

1. Basic Properties of Stars



This is the Sun during a total eclipse. The Sun, our closest star, is very much representative of the objects that we will study during this module, namely stars. Much of the knowledge that we currently have about stars has been derived by studying the Sun. In this module, we start by introducing a number of important observations that we have of stars, and that are important to understand their structure. We then continue by dedicating the largest part of the module to studying the interiors of stars. We end the module by talking about stellar structure: first in a schematic way, and then more phenomenological.

1.1 Basic Properties of Stars

- a) Luminosity L
- b) Effective Temperature T_{eff}
- c) Chemical Composition X, Y, Z
- d) Radius

e) Mass

f) Age

With these properties of stars we should be able to describe most of their formation, evolution and death. In the next part of this chapter, we will define these quantities, explain how they are being obtained, and give representative values for the only star in our Solar System: the Sun.

For the Sun we have:

$$L_{\text{TM}} = 4 \cdot 10^{26} \text{ W}$$

$$M_{\text{TM}} = 2 \cdot 10^{30} \text{ kg}$$

$$R_{\text{TM}} = 7 \cdot 10^5 \text{ km}$$

$$(T \sim 6000\text{K})$$

$$(\text{Age} \sim 10^{10} \text{ y})$$

For other stars we have the following range:

$$10^{-4} \leq \frac{L}{L_{\text{Sun}}} \leq 10^6$$

$$\frac{1}{3} \leq \frac{T_{\text{eff}}}{T_{\text{eff}, \text{Sun}}} \leq 3$$

$$10^{-2} \leq \frac{R}{R_{\text{Sun}}} \leq 10^3$$

$$10^{-1} \leq \frac{M}{M_{\text{Sun}}} \leq 10^2$$

1.2 Measured Quantities

a) Luminosity

We cannot directly measure the luminosity of a star, since the amount of light detected, the flux, depends on the distance.

$$\text{Energy Flux} = \frac{\text{Power output}}{\text{Area}} = \frac{L}{4 \pi d^2} = b$$

Here d is the distance of the stars.

For example, on a photograph one can see many thousands of stars, some brighter than others. If a star is brighter than another on such a plate, it doesn't mean that its luminosity is larger. It can just as well mean that it is nearer.



The quantity $b = \frac{L}{4\pi d^2}$ is sometimes called the bolometric flux, i.e., the flux measured by a detector with perfect sensitivity at all wavelengths.

We define:

$$m_{bol} \equiv -2.5 \log_{10} \left(\frac{L}{4\pi d^2} \right) + cst.$$

m_{bol} is here called the apparent bolometric magnitude.

The constant in the past was determined by requiring that the stars Vega (α Lyr) had a bolometric magnitude 0.00. Later this has been changed by defining this zero-point based on a number of nearby stars.

This definition is equivalent with:

$$m_{bol} = -2.5 \log(L) + 5 \log(d) + cst.$$

(From now on we use \log for \log_{10})

The definition implies that a brighter object has a lower magnitude. Also, an object that is a factor 10 further away is 5 magnitudes fainter. As an example, $m_{bol, Sun} = -27.1$

Exercise: Assume that the moon reflects all the light from the Sun, what is then its apparent magnitude?

Answer: The brightness of the Sun is: $b_{Sun} = \frac{L_{Sun}}{4 \pi d_{Sun}^2}$

This means that $b_{moon} = b_{Sun} \frac{\pi R_{moon}^2}{4 \pi d_{moon}^2}$, where R_{moon} = the radius of the moon, and

d_{moon} is the distance from the Earth to the Moon.

We have then:

$$\frac{b_{moon}}{b_{Sun}} = \frac{R_{moon}^2}{4 \pi d_{moon}^2} \sim \frac{(10^3 \text{ km})^2}{4 \pi (3.8 \cdot 10^5)^2} \sim 10^{-5}$$

Now convert this brightness ratio into a magnitude difference:

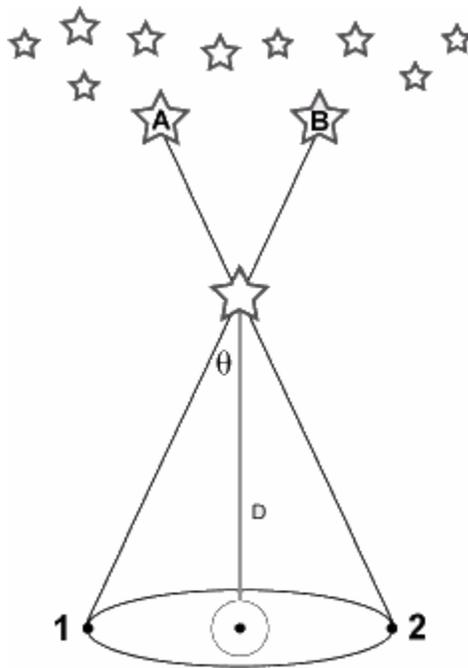
$$m_{moon} - m_{Sun} = -2.5 \log\left(\frac{b_{Moon}}{b_{Sun}}\right) = 12.5$$

Since $m_{bol, Sun} = -27.1$ we have $m_{bol, moon} \sim -14.6$

To explain why we use these *magnitudes* one has to go back to the Greeks, who classified stars in 6 brightness classes 1 (brightest) to 6 (faintest). Since they had been defined by eye, and the eye has a logarithmic response, it turned out that a magnitude corresponds to about a factor 2.5 in brightness. The new system has been defined in such a way as to be able to more or less produce the old star catalogues.

To be able to determine luminosities we need to know the distance of stars. This is a complicated problem. More details about how distances can be measured can be found here. For nearby stars an accurate method for the determination of distances is the *parallax method*.

The most direct method to measure the distance to nearby stars is through the use of parallax. The Earth's motion around the Sun every year produces a very small shift in nearby star's position in the sky compared to distant, background stars. This shift is



Parallax

always less than one arcsec for any star, which is very small (where a circle is 360 degrees, one degree is 60 arcminutes, one arcminute is 60 arcsecs).

The parallax of a star is used to calculate its distance. The angle of the parallax, θ , in arcsecs is equal to the distance, D by the following formula:

$$D = \frac{1}{\theta}$$

here the distance, D is measured in parsecs, where one parsec equals 3.26 light-years (the distance light travels in one year).

An object for which the parallax is 1 arcsec is by definition at the distance of 1 pc. This corresponds to 206265 AU or $3.086 \cdot 10^{16}$ m. Based on the luminosity, we define the distance-independent *absolute magnitude* in the following way:

$$M_{bol} \equiv -2.5 \log\left(\frac{L}{10 pc}\right) + cst$$

, where the constant is determined by requiring that the star

Vega (α Lyr) has absolute bolometric magnitude 0.

Since we also had:

$$m_{bol} = -2.5 \log\left(\frac{L}{4\pi d^2}\right) + cst$$

we have

$$m_{bol} - M_{bol} = -5 + 5 \log d$$

This is an important equation, which makes it easy to obtain absolute magnitudes from apparent magnitudes.

Question: What is the absolute magnitude of the Sun?

Answer: We had that $m_{bol} = -27.1$

$$m_{bol} - M_{bol} = -5 + 5 \log d = 5 \log \left(\frac{d}{10 \text{ pc}} \right) = 5 \log (3 \cdot 10^{-7})$$

So $M_{bol} = +5.5$

b. Effective Temperature

Stars are radiating objects, and a first approximation to their spectra can be a black-body curve:

$$B_\nu(T) = \frac{2h\nu^3}{c^2} \frac{1}{e^{\frac{h\nu}{kT}} - 1}, \text{ or } b_\lambda(T) = \frac{2hc^2}{\lambda^5} \frac{1}{e^{\frac{hc}{\lambda kT}} - 1}$$

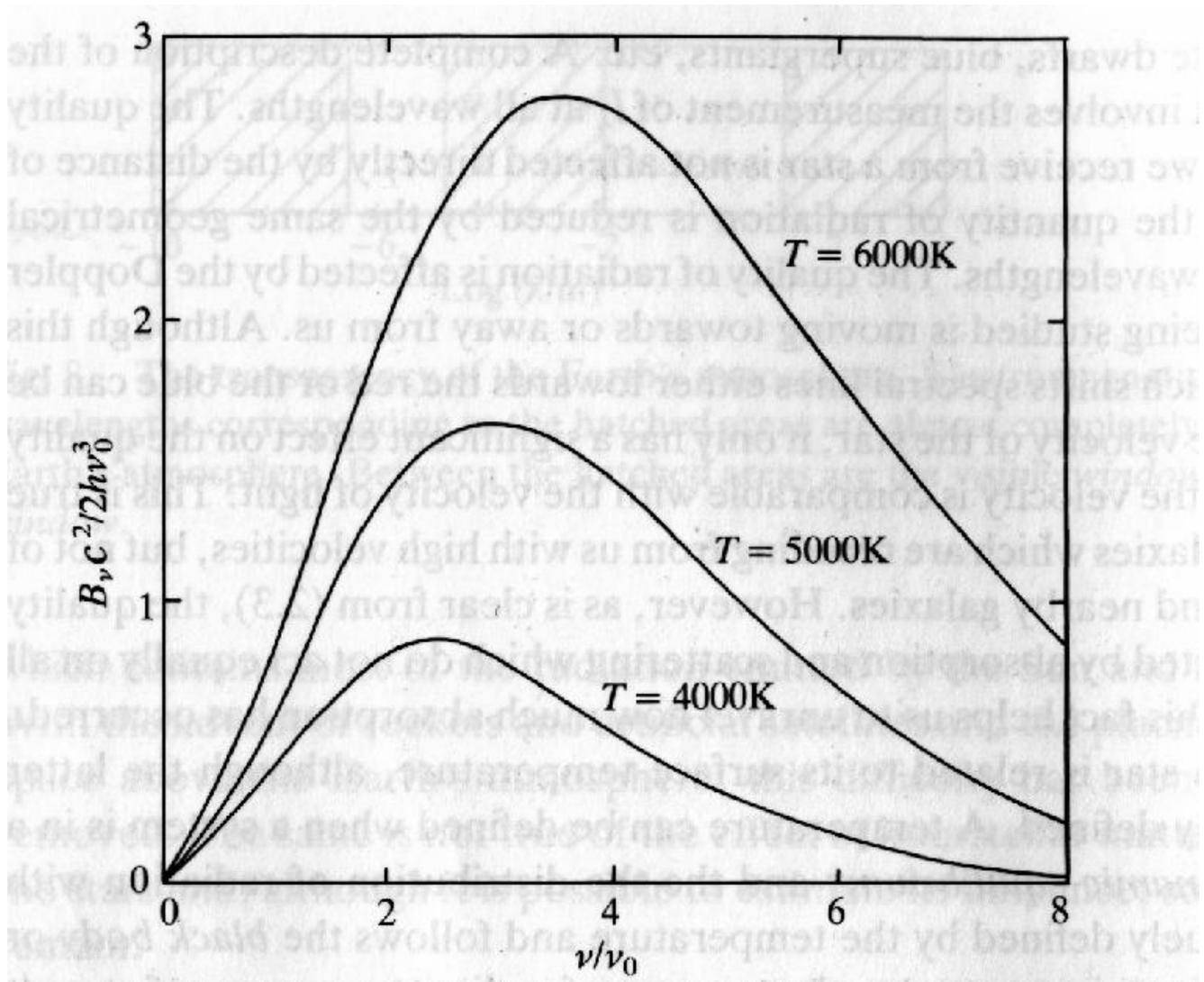
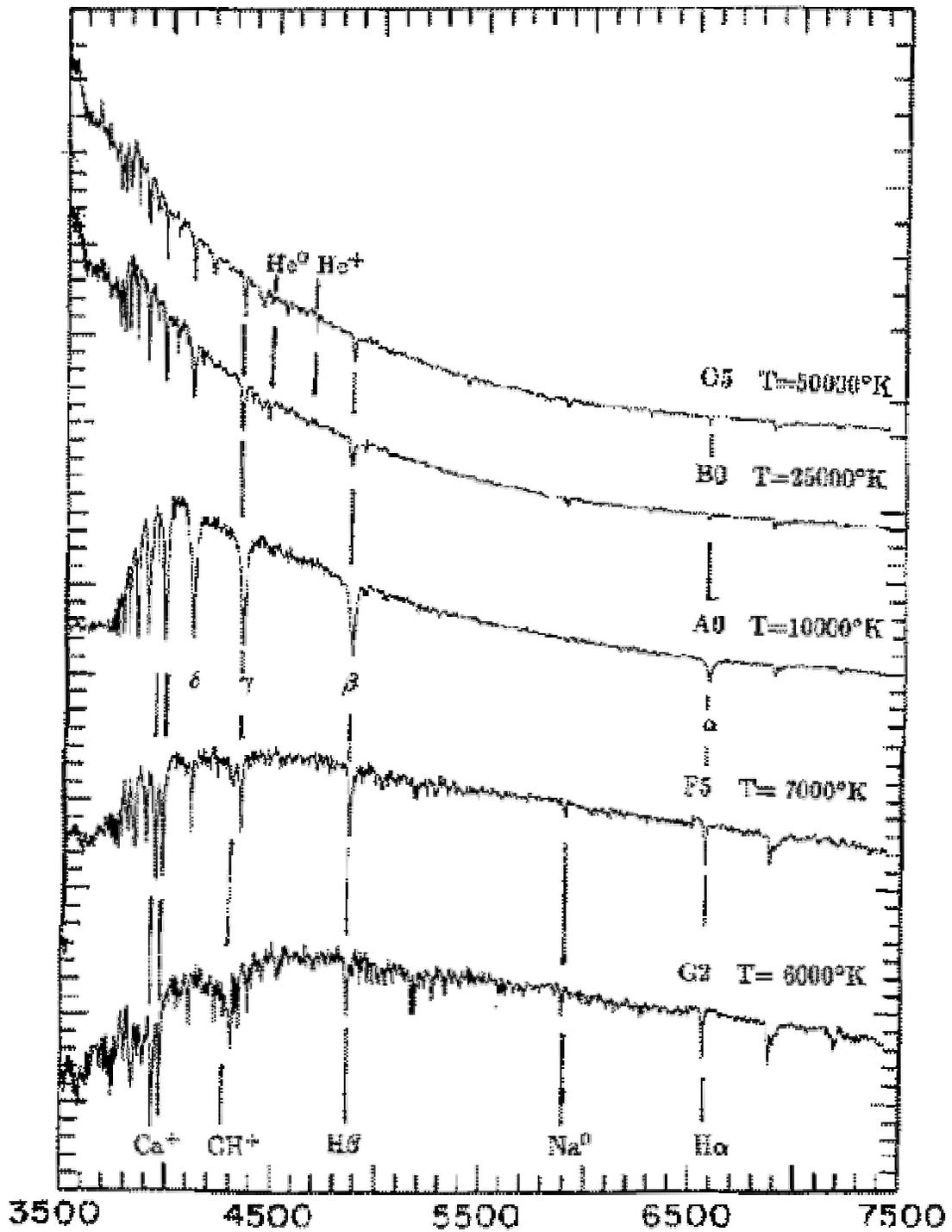


Fig. 1: Some blackbody curves.

In fact, for many stars the blackbody approximation is not such a good one, since their spectrum contain absorption and emission lines, and might have several thermal components in their spectrum (e.g, photosphere, corona, etc.).

Here some spectra of stars:



We now define the *effective temperature* of a star as the black body temperature of an object with the same radius that radiates the same amount of energy.
 So, for a blackbody with radius R and luminosity L the temperature would be defined by

$$L = 4 \pi R^2 \sigma T^4$$

For a star with the same radius and luminosity we define the effective temperature as

$$L = 4 \pi R^2 \sigma T_{eff}^4$$

This is Stefan-Boltzmann's law. Here σ is the constant of Stefan-Boltzmann, which is about $5.6710^{-8} \text{ W m}^{-2} \text{ K}^{-4}$

Since stars are often very faint, one cannot easily measure monochromatic fluxes. For that reason one often uses *colours* in astronomy.

c) Colours

Colours are defined by comparing power outputs over different parts of the spectrum. Let the bolometric luminosity be

$$L_{bol} = \int_0^{\infty} L(\lambda) d\lambda$$

In practise one always measure the flux in a certain bandpass. So for a filter with bandpass $f(\lambda)$ we have

$$L_f = \int_0^{\infty} L(\lambda) f(\lambda) d\lambda$$

Astronomers have defined several filter systems consisting of a number of filters, that can be used to quickly measure observational parameters of stars. For example, the most commonly used filter system, Johnson (1966)'s UBV system contains 3 filters:

Filter	Central wavelength	Filter Width
U	360 nm	70 nm
B	440 nm	100 nm
V	550 nm	90 nm

By subtracting 2 bands one obtains the slope of the spectrum between two characteristic wavelengths. This slope correlates with temperature, so to first order a *colour* (e.g. B-V or U-B) is a good measure of the temperature of a star. A second colour then gives information about other parameters, like chemical abundance.

The U-band flux is defined by

$$U = -2.5 \log \left(\int_0^{\infty} \frac{f_U(\lambda) L(\lambda) d\lambda}{4 \pi d^2} \right) + cst_U$$

for a different band, say V, we have

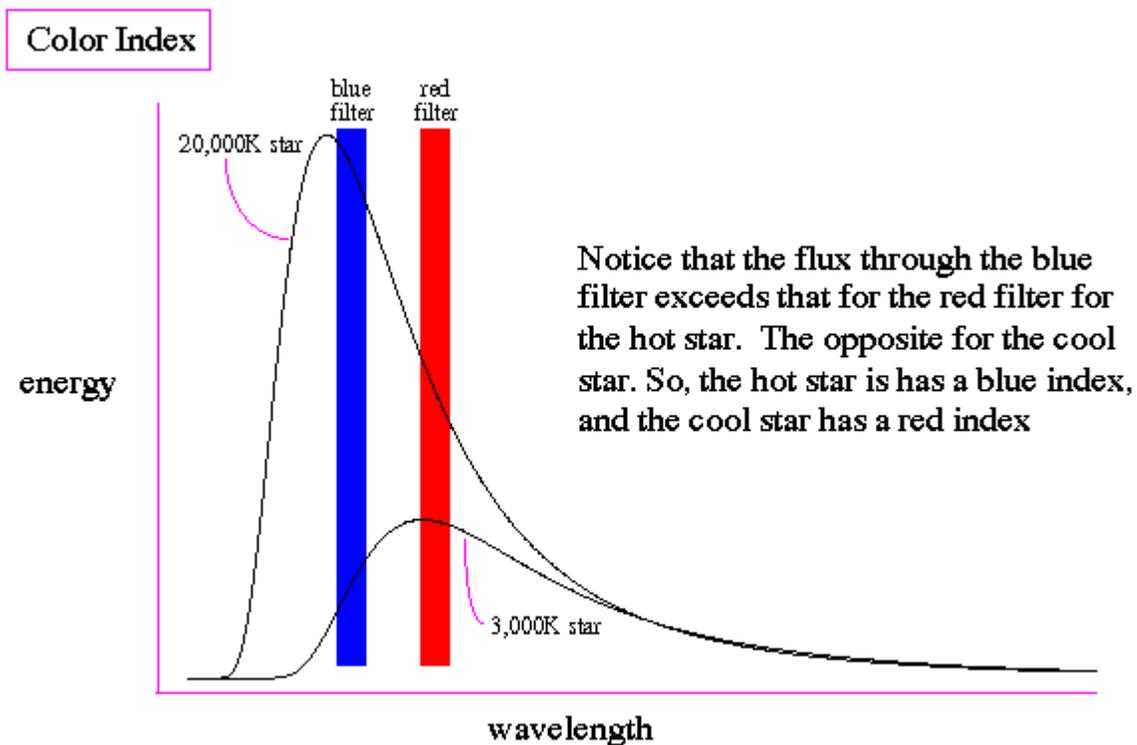
$$V = -2.5 \log \left(\int_0^\infty \frac{f_V(\lambda) L(\lambda) d\lambda}{4\pi d^2} \right) + cst_V$$

so $U - V = cst - 2.5 \log \left(\frac{\int_0^\infty L(\lambda) f_U(\lambda) d\lambda}{\int_0^\infty L(\lambda) f_V(\lambda) d\lambda} \right)$ is independent of distance d.

Roughly, we can approximate B-V by:

$$B - V = \frac{7000 K}{T_{eff}} - 0.56$$

(showing that cooler, redder stars have a larger B-V).



One often tries to determine the luminosity of a star by measuring the magnitude in a certain band, most often V. The error that one makes in estimating the total bolometric magnitude in this way depends on the distance of the effective wavelength of the star from the centre of the V-band. One defines the bolometric correction B.C as

$$B.C. \equiv M_{bol} - M_V$$

Bolometric corrections are often tabulated for all kinds of stars, making it possible to easily determine bolometric magnitudes.

d) The Electromagnetic Spectrum

The spectrum of a star is basically the spectrum of a blackbody with a temperature between 2000 and 20000 K, with absorption lines due to transitions near the surface of the star (this zone is called the photosphere) as well as some emission lines due to transitions of ions above the photosphere. Since hydrogen is by far the most abundant element, transitions between different states of hydrogen are very important (see later).

For hydrogen, with one proton and one electron, we have the following energy levels (see the first year's lectures):

$$E_n = \frac{\mu e^4}{2\hbar^2} \frac{1}{n^2} = -\frac{13.6}{n^2} eV$$

Here μ is the mass of the electron, and e its charge.

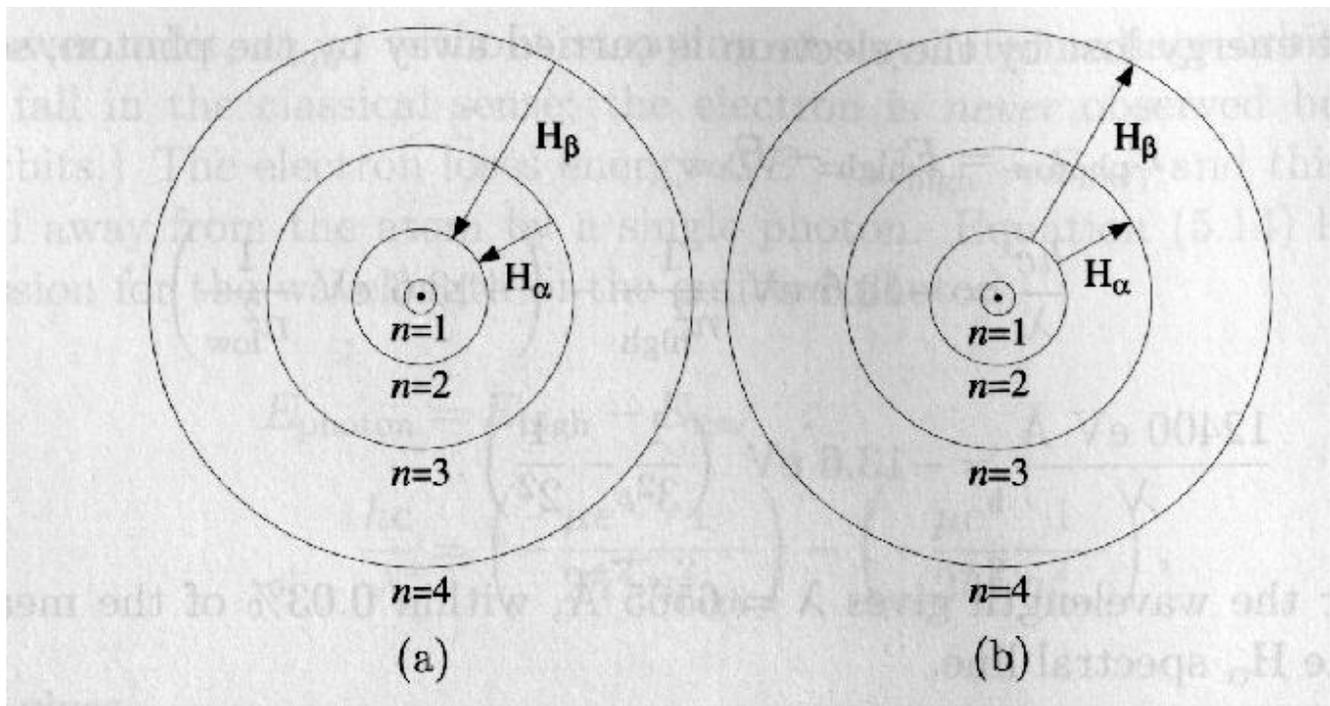


Figure: Balmer absorption lines produced by the Bohr atom. (a) absorption lines, and (b) emission lines.

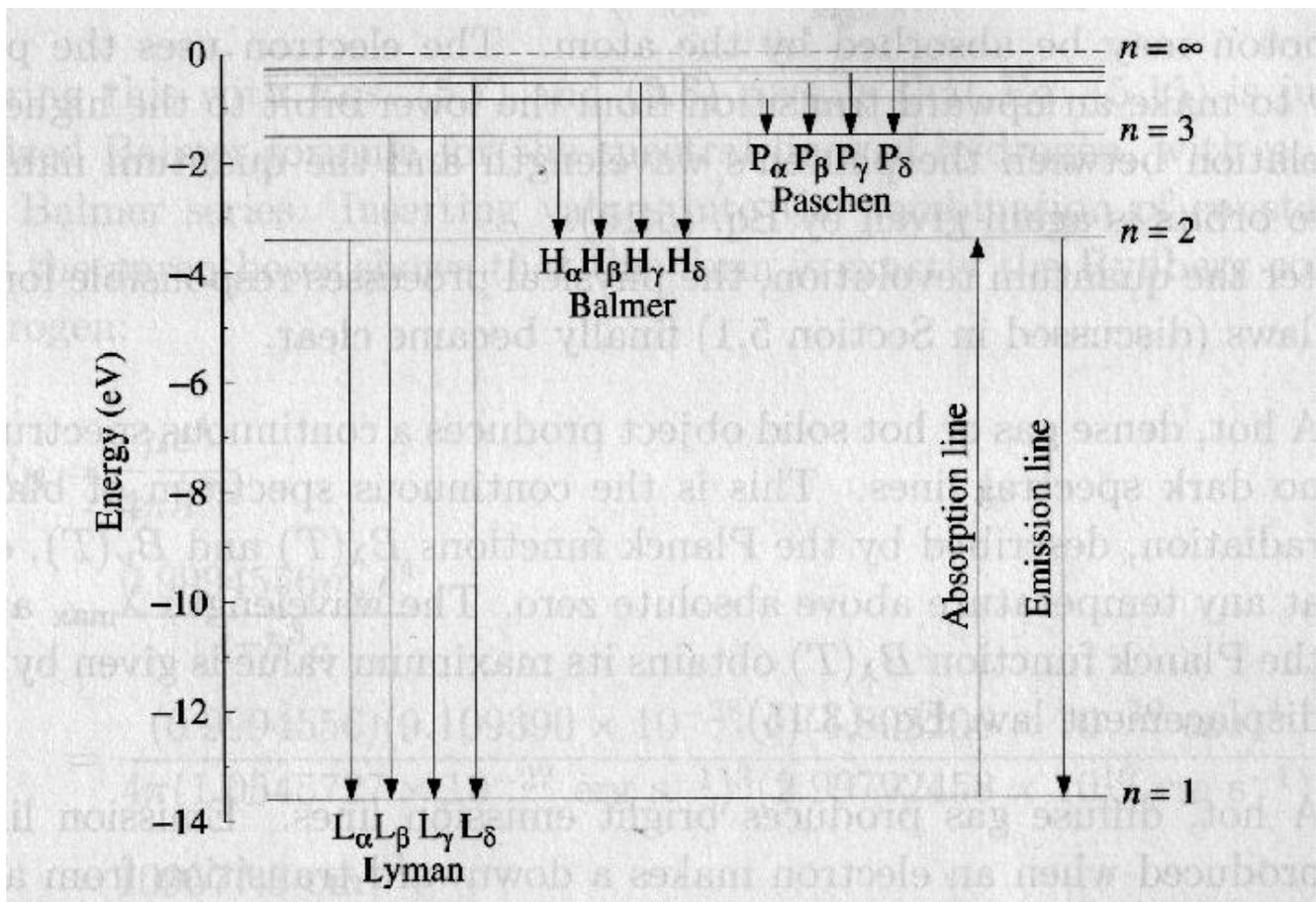


Figure: Energy level diagram for the Balmer lines.

The transitions from E_n to E_1 are called the *Lyman* series. The transitions to E_2 are called the *Balmer* series, while the transitions to E_3 are called the *Paschen* series. For example, the transition from E_3 to E_2 is called $H\alpha$, and is the first line of the Balmer series.

e) Stellar Masses

Since stars are so far away, it is impossible to measure their masses directly. Instead, we look for binary star systems and use Newton's law of gravity to measure their masses.

Two stars in a binary system are bound by gravity and revolve around a common center of mass. Kepler's 3rd law of planetary motion can be used to determine the sum of the mass of the binary stars if the distance between each other and their orbital period is known.

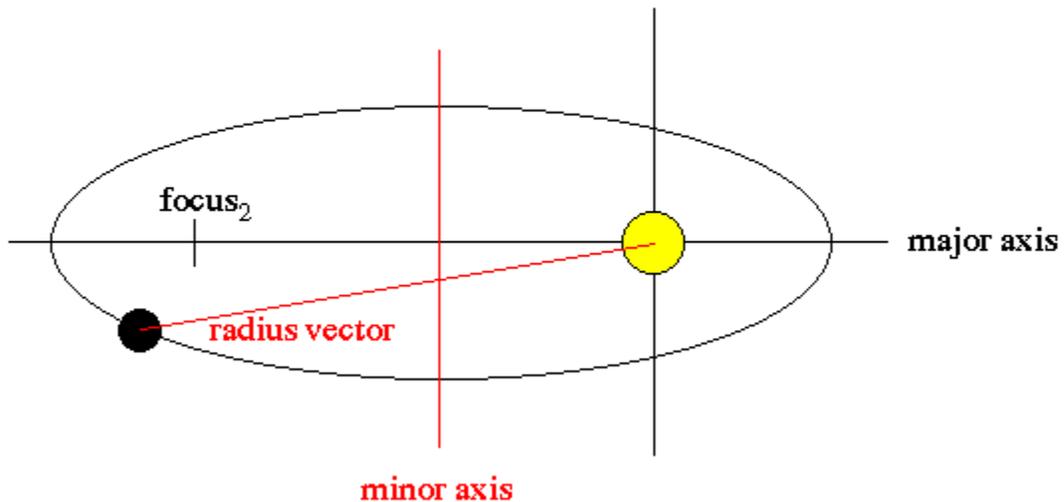
If two stars orbit each other at large separations, they evolve independently and are called a wide pair. If the two stars are close enough to transfer matter by tidal forces,

then they are called a close or contact pair.

Binary stars obey Kepler's Laws of Planetary Motion, of which there are three.

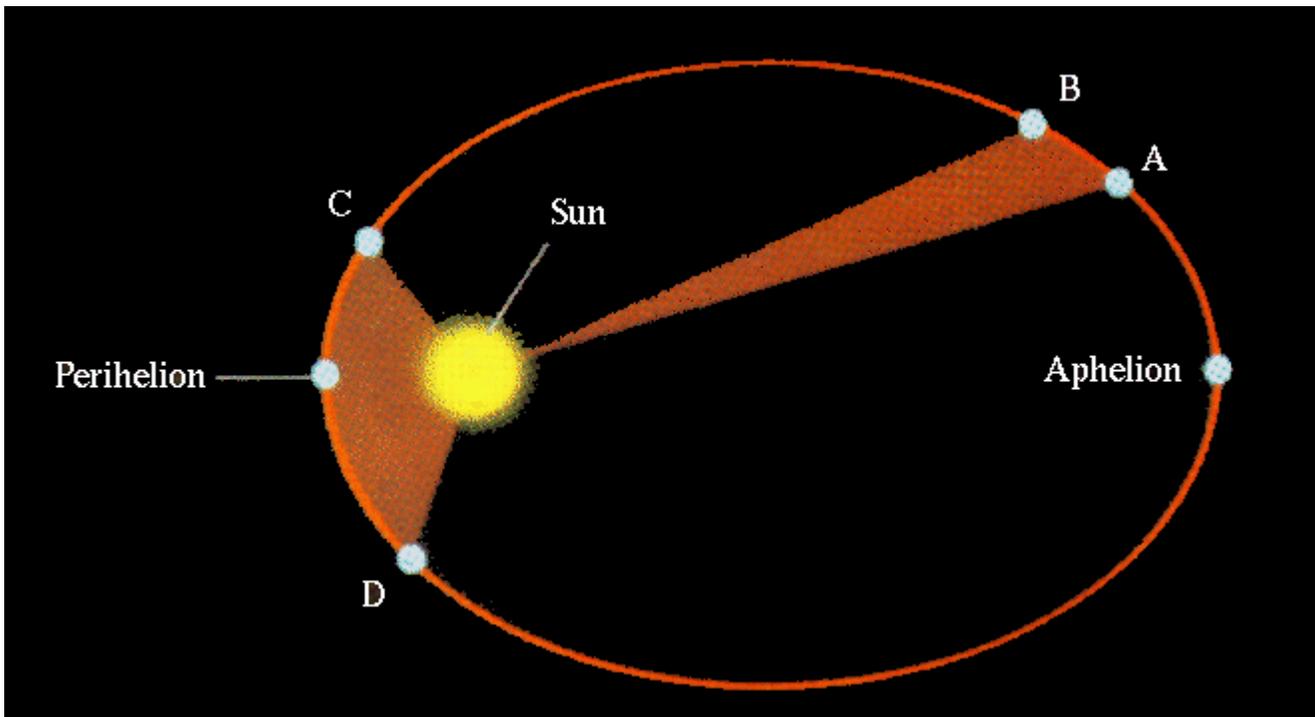
1st law (law of elliptic orbits): Each star or planet moves in an elliptical orbit with the center of mass at one focus.

Ellipse = member of a family of mathematical curves called conic sections



Ellipses that are highly flattened are called highly eccentric. Ellipses that are close to a circle have low eccentricity.

2nd law (law of equal areas): a line between one star and the other (called the radius vector) sweeps out equal areas in equal times



This law means that objects travel fastest at the low point of their orbits, and travel slowest at the high point of their orbits.

3rd law (law of harmonics): The square of a star or planet's orbital period is proportional to its mean distance from the center of mass cubed

It is this last law that allows us to determine the mass of the binary star system (note only the sum of the two masses):

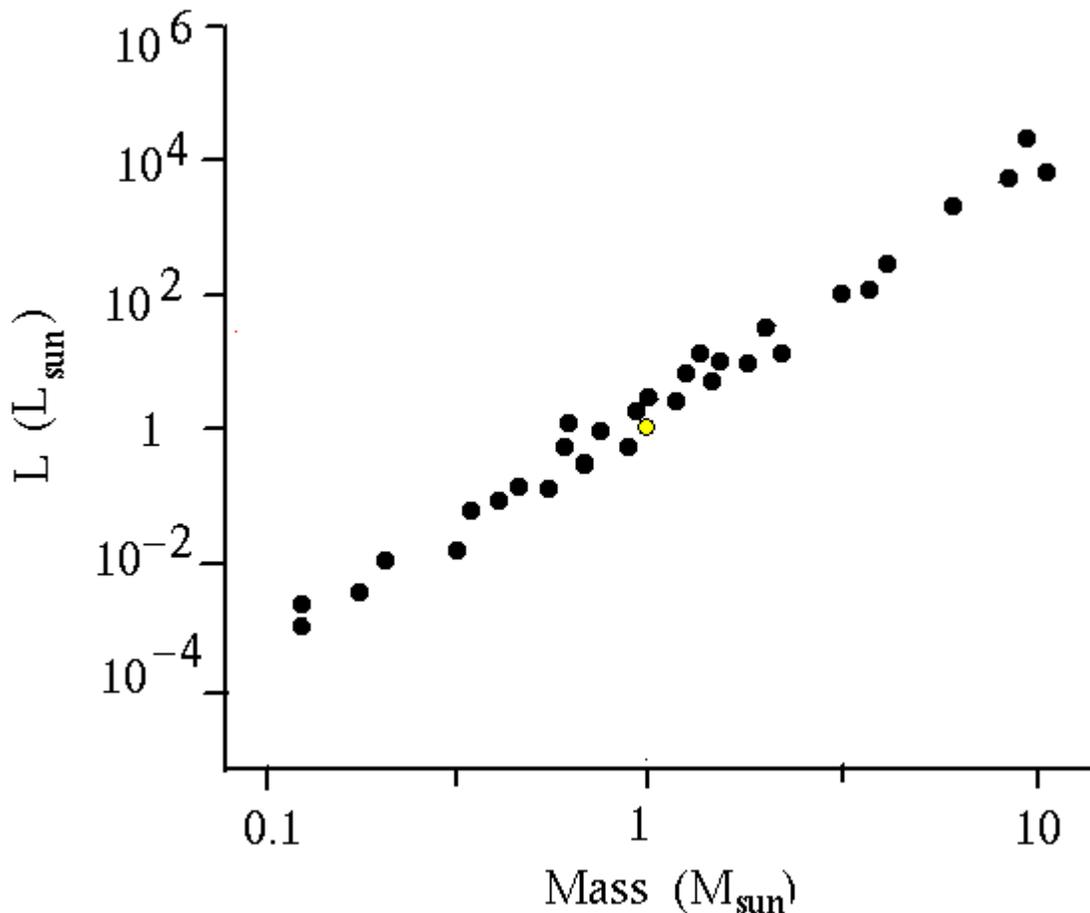
Kepler's 3rd law states the square of a planet or stars orbital period is proportional to its mean distance from each other such that:

$$r^3 = kP^2$$

where P is the orbital period in years and r is the distance between each other in Astronomical Units (the distance from the Earth to the Sun). The constant, k , is derived from Newton's law of gravity to be the sum of the masses of the stars, $M_1 + M_2$, in units of solar mass. So the full equation becomes:

$$M_1 + M_2 = \frac{r^3}{P^2}$$

When you plot the mass of a star versus its absolute luminosity, one finds a correlation between the two quantities shown below.



This relationship is called the [mass-luminosity relation](#) for stars, and it indicates that the mass of a star controls the rate of energy production, which is thermonuclear fusion in the star's core. The rate of energy generation, in turn, uniquely determines the star's total luminosity. Note that this relation only applies to stars before they evolve into giant stars (those stars which burn hydrogen in their core).

The above diagram is for stars near the Sun, a small sample. When we examine all the stars in our Galaxy we find that they range in mass from about 0.08 to 100 times the mass of the Sun. The lower mass limit is set by the internal pressures and temperatures needed to start thermonuclear fusion (protostars too low in mass never beginning fusion and do not become stars). The upper limit is set by the fact that stars of mass higher than 100 solar masses become unstable and explode. Notice also that these range of masses corresponds to a luminosity range from 0.0001 to 10^5 solar luminosities.

f) Stellar Radii

Stellar radii are small. Their diameter cannot be determined from direct imaging observations. For example, a star like the Sun at the distance of the brightest star, Proxima Centauri, would only have an angular size of $6 \cdot 10^{-3}$ arcsec.

There are 3 often used methods to determine stellar radii:

1. Using Stefan-Boltzmann's law.

Stefan-Boltzmann's law relates the energy emitted by an object to its temperature. However, this is for point sources, an idealized case. Real objects have size, which means surface area. Large objects cool faster (emit energy faster) than small objects. For stars, the luminosity, temperature and radius are linked by an expansion of Stefan-Boltzmann's law:

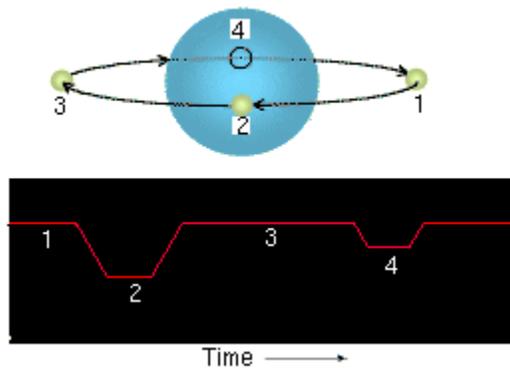
$$L = 4\pi R^2 \sigma T^4$$

This equation can be expressed in terms of solar units such that:

$$\frac{L}{L_{\odot}} = \left[\frac{R}{R_{\odot}} \right]^2 \left[\frac{T}{T_{\odot}} \right]^4$$

This needs knowledge of d , to get the Luminosity, and the electromagnetic spectrum.

2. Using eclipsing binaries.



For an eclipsing binary, one gets the radius of the stars by measuring the time that the

star is being eclipsed by the other. To do this, we also need the distance.

3. Using interferometric methods

One uses an interferometer (i.e. 2 radio telescopes close together) to measure the radius of a star.

g) Chemical Composition and Spectral Types

From absorption and emission lines the chemical composition in the photosphere of the star can be determined. We define

X = the relative mass fraction of hydrogen

Y = the relative mass fraction of helium

Z = the relative mass fraction of all other elements.

For the Sun $X=0.75$, $Y=0.23$ and $Z=0.02$

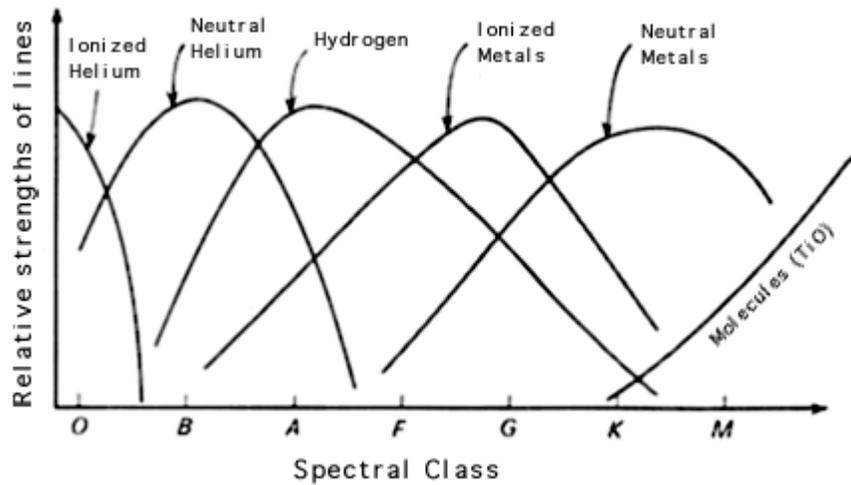
Spectral Types

Stars are divided into a series of spectral types based on the appearance of their absorption spectra. Some stars have a strong signature of hydrogen (O and B stars), others have weak hydrogen lines, but strong lines of calcium and magnesium (G and K stars). After years of cataloging stars, they were divided into 7 basic classes: O, B, A, F, G, K and M. Note that the spectra classes are also divisions of temperature such that O stars are hot, M stars are cool.

Between the classes there were 10 subdivisions numbered 0 to 9. For example, our Sun is a G2 star. Sirius, a hot blue star, is type B3.

Why do some stars have strong lines of hydrogen, others strong lines of calcium? The answer was not composition (all stars are 95% hydrogen) but rather surface temperature. As temperature increases, electrons are kicked up to higher levels (remember the Bohr model) by collisions with other atoms. Large atoms have more kinetic energy, and their electrons are excited first, followed by lower mass atoms.

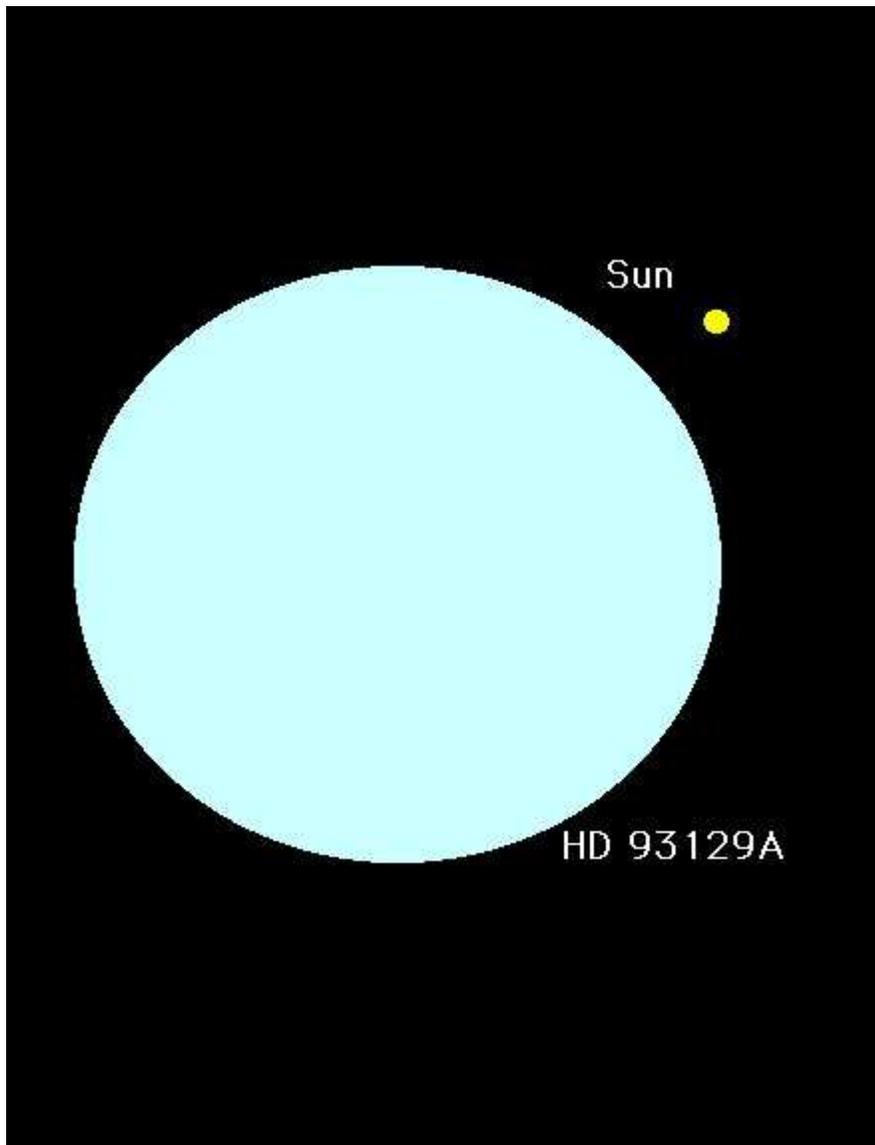
If the collision is strong enough (high temperatures) then the electron is knocked off the atom and we say the atom is ionized. So as we go from low temperatures in stars (couple 1,000K) we see heavy atoms, like calcium and magnesium, in the stars spectrum. As the temperature increases, we see lighter atoms, such as hydrogen (the heavier atoms are all ionized by this point and have no electrons to produce absorption lines).



As we will see later, hotter stars are also more massive stars (more energy burned in the core). So the spectral classes of stars is actually a range of masses, temperatures, sizes and luminosity. For normal stars (called main sequence stars) the following table gives their properties:

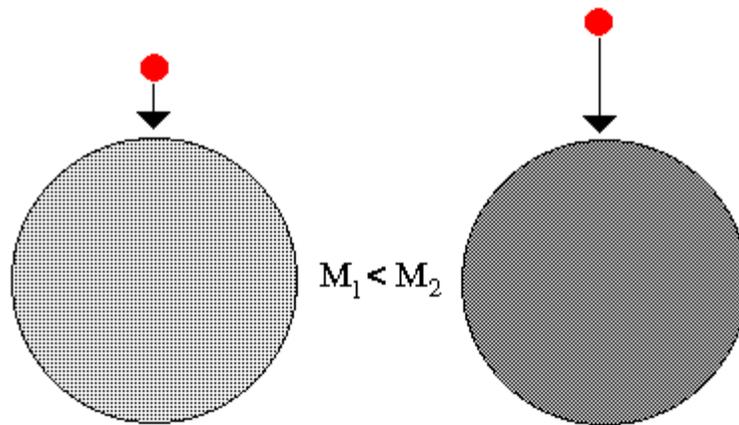
type	Mass	Temp	Radius	Lum (Sun=1)
O	60.0	50,000	15.0	1,400,000
B	18.0	28,000	7.0	20,000
A	3.2	10,000	2.5	80
F	1.7	7,400	1.3	6
G	1.1	6,000	1.1	1.2
K	0.8	4,900	0.9	0.4
M	0.3	3,000	0.4	0.04

So our Sun is a fairly middle-of-the-road G2 star: A B star is much larger, brighter and hotter. An example is HD93129A shown below:

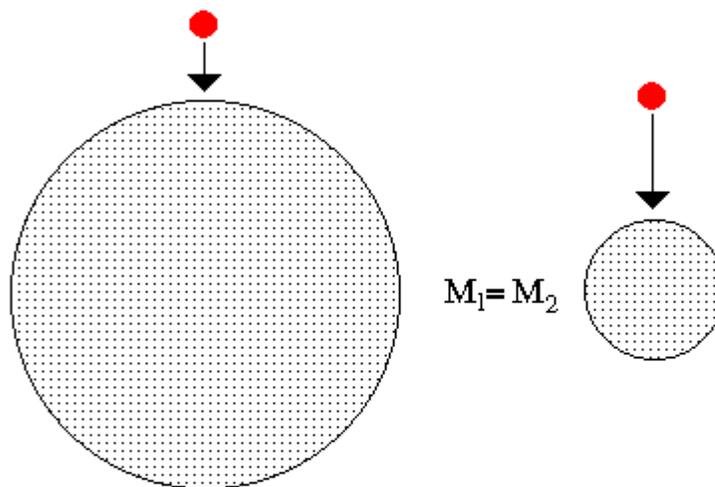


Luminosity Classes:

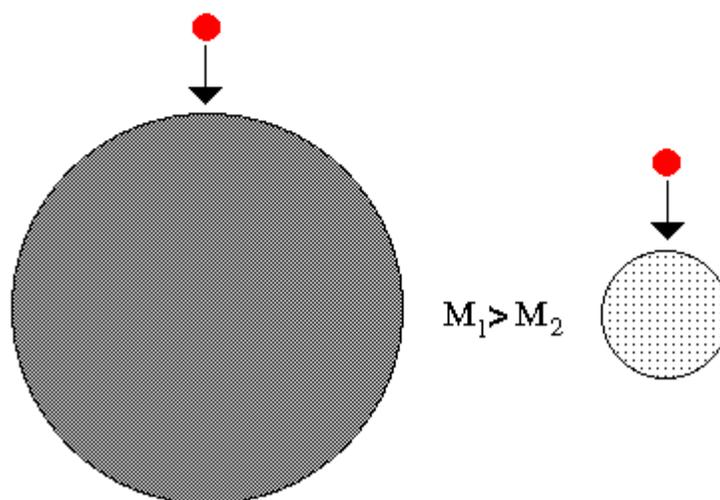
Closer examination of the spectra of stars shows that there are small changes in the patterns of the atoms that indicate that stars can be separated by size called luminosity classes. The strength of a spectra line is determined by what percentage of that element is ionized. An atom that is ionized has had all its electrons stripped off and can produce no absorption of photons. At low densities, collisions between atoms are rare and they are not ionized. At higher densities, more and more of the atoms of a particular element become ionized, and the spectral lines become weak. One way to increase density at the surface of a star is by increasing surface gravity. The strength of gravity at the surface of a star is determined by its mass and its radius (remember escape velocity). For two stars of the same mass, but different sizes, the larger star has a lower surface gravity = lower density = less ionization = stronger spectral lines.



as the mass of the star increases, the surface gravity increases



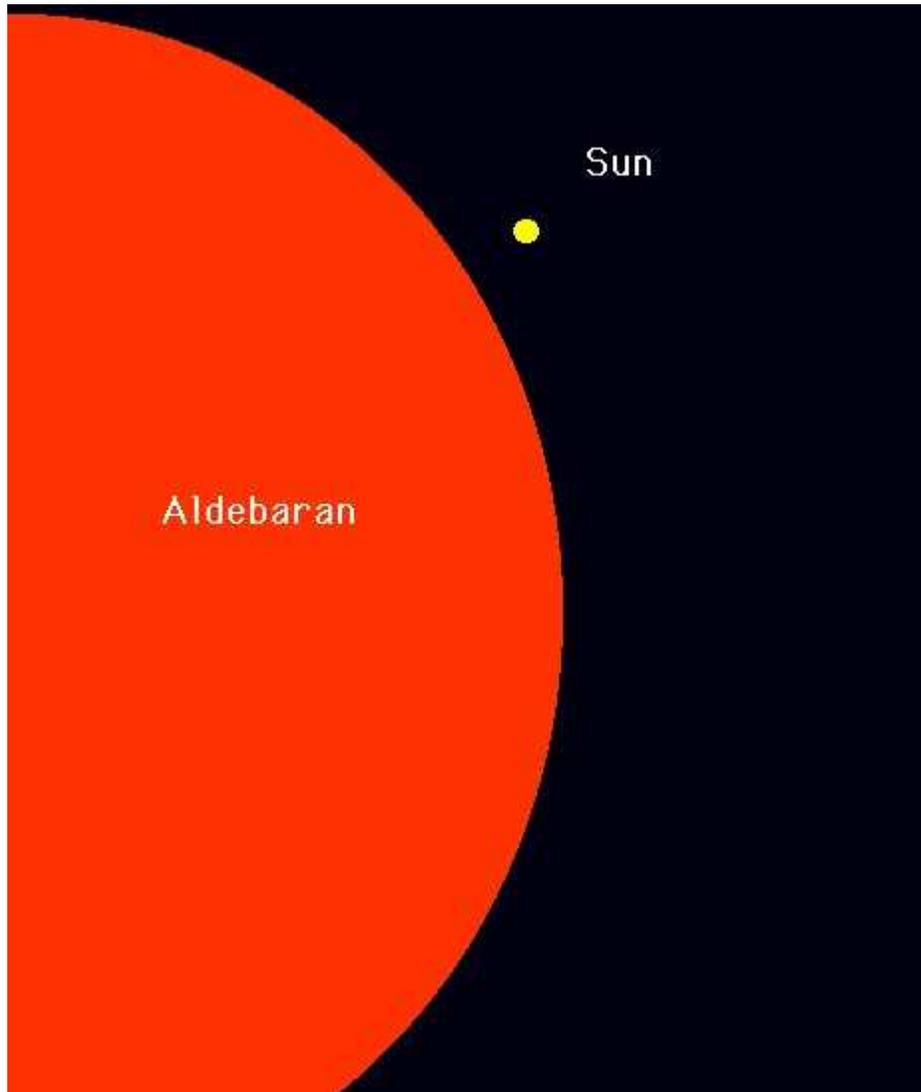
as the radius of the star decreases, the surface gravity increases



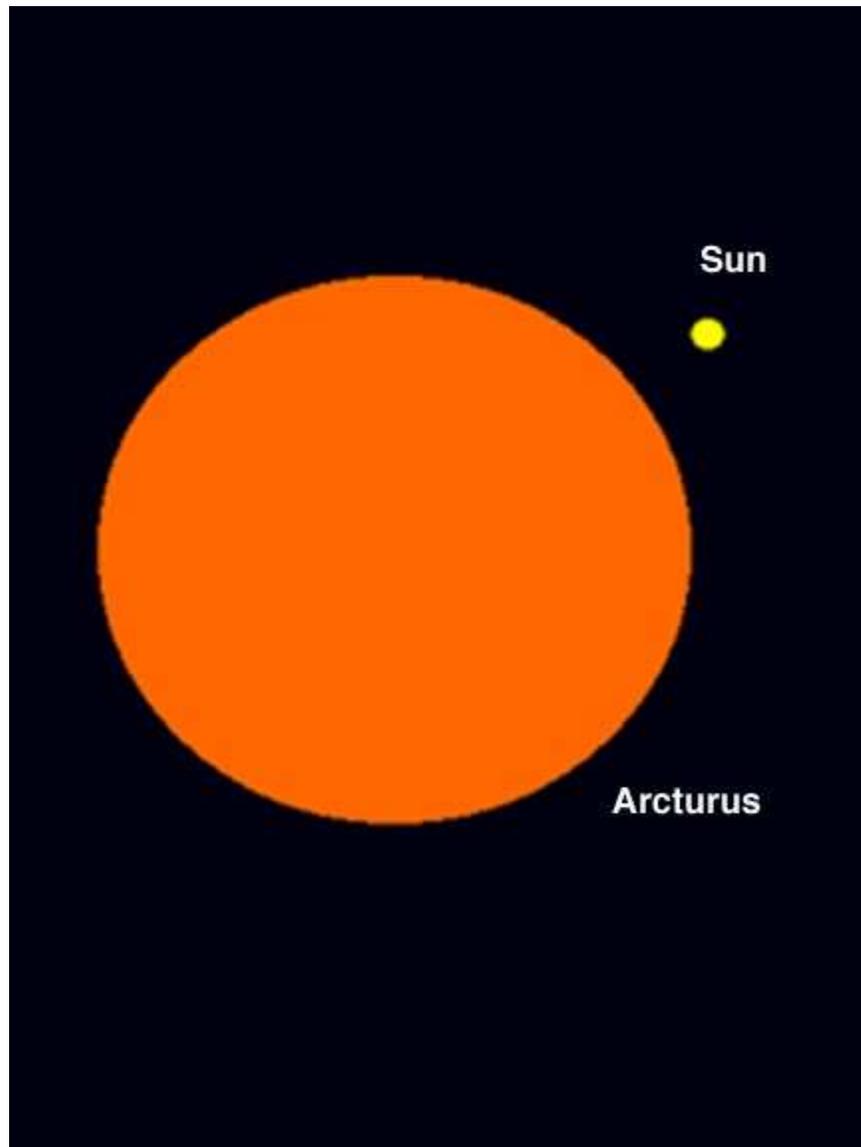
the surface gravity of a large, massive star can equal the surface gravity of the small, less massive star.

This was applied to all stars and it was found that stars divide into five luminosity classes: I, II, III, IV and V. Stars of type I and II are called supergiants, being very large (low surface gravity), stars of type III and IV are called giant stars. Stars of type V are called dwarfs. The Sun is a G2 V type stars.

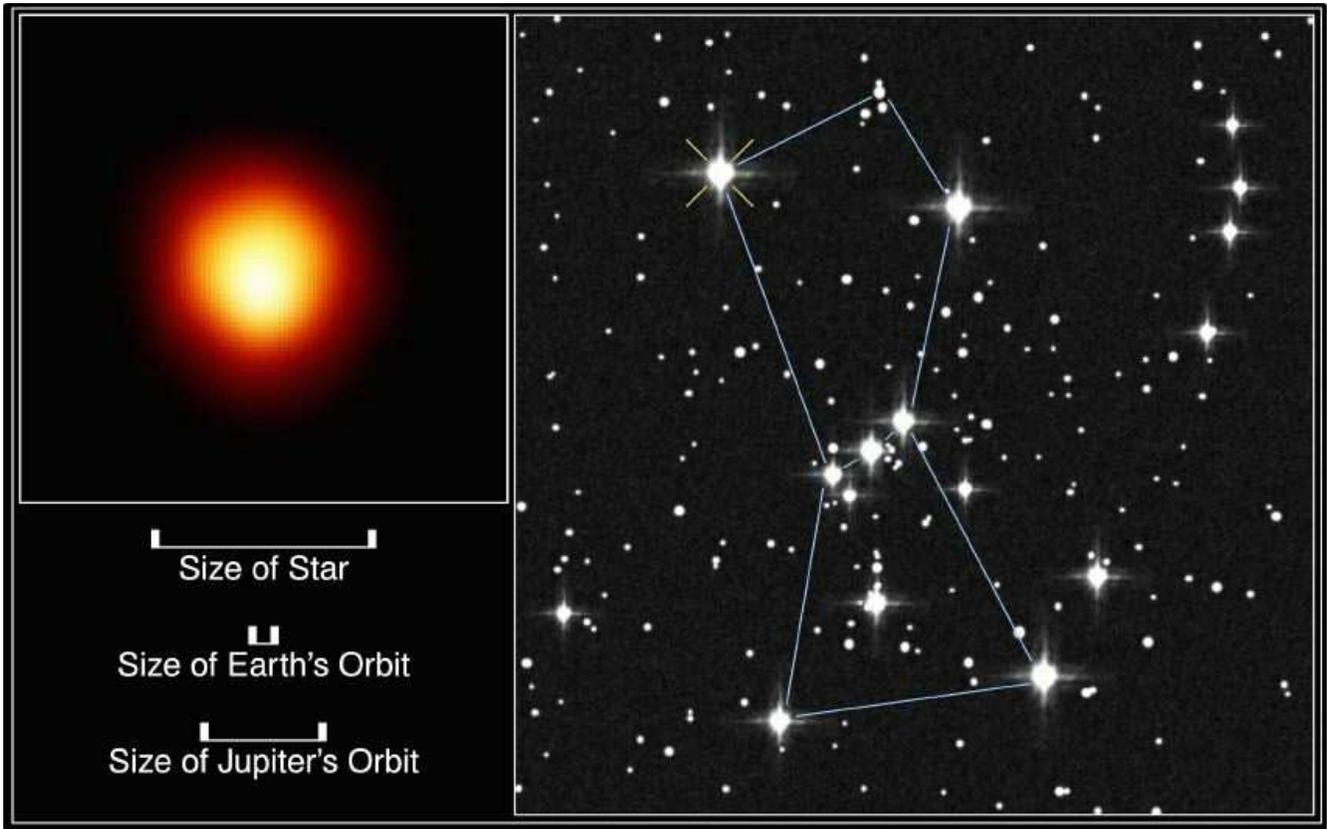
So now we have a range of stellar colors and sizes. For example, Aldebaran is a red supergiant star:



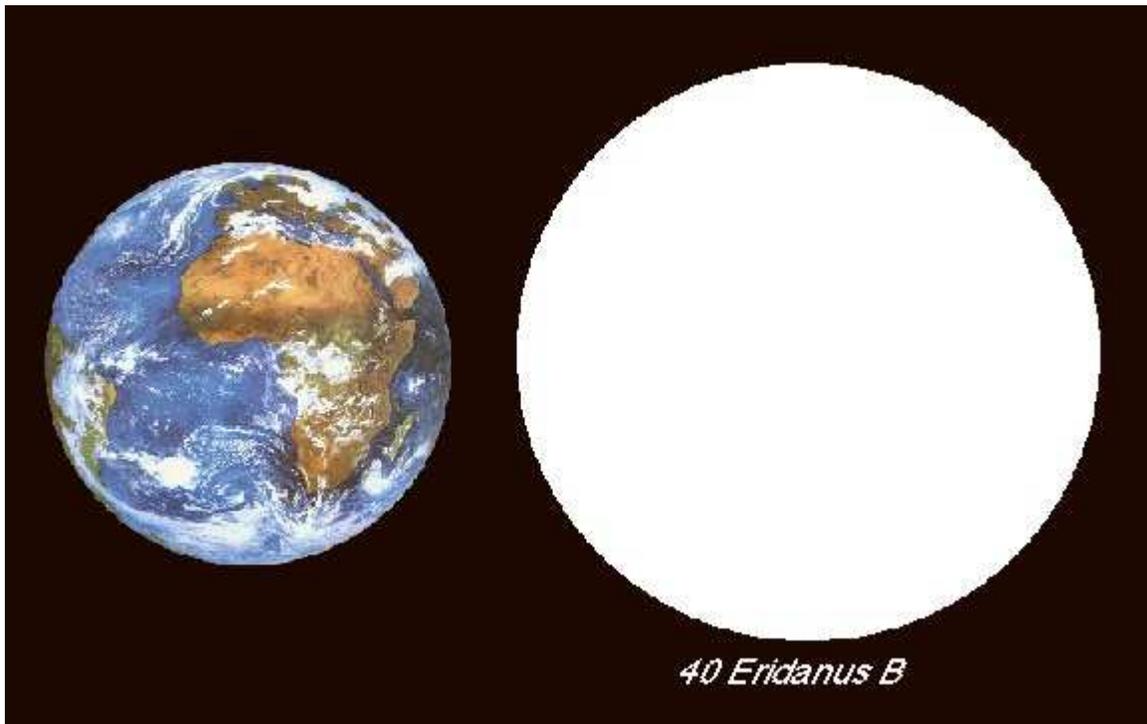
Arcturus is an orange giant star:



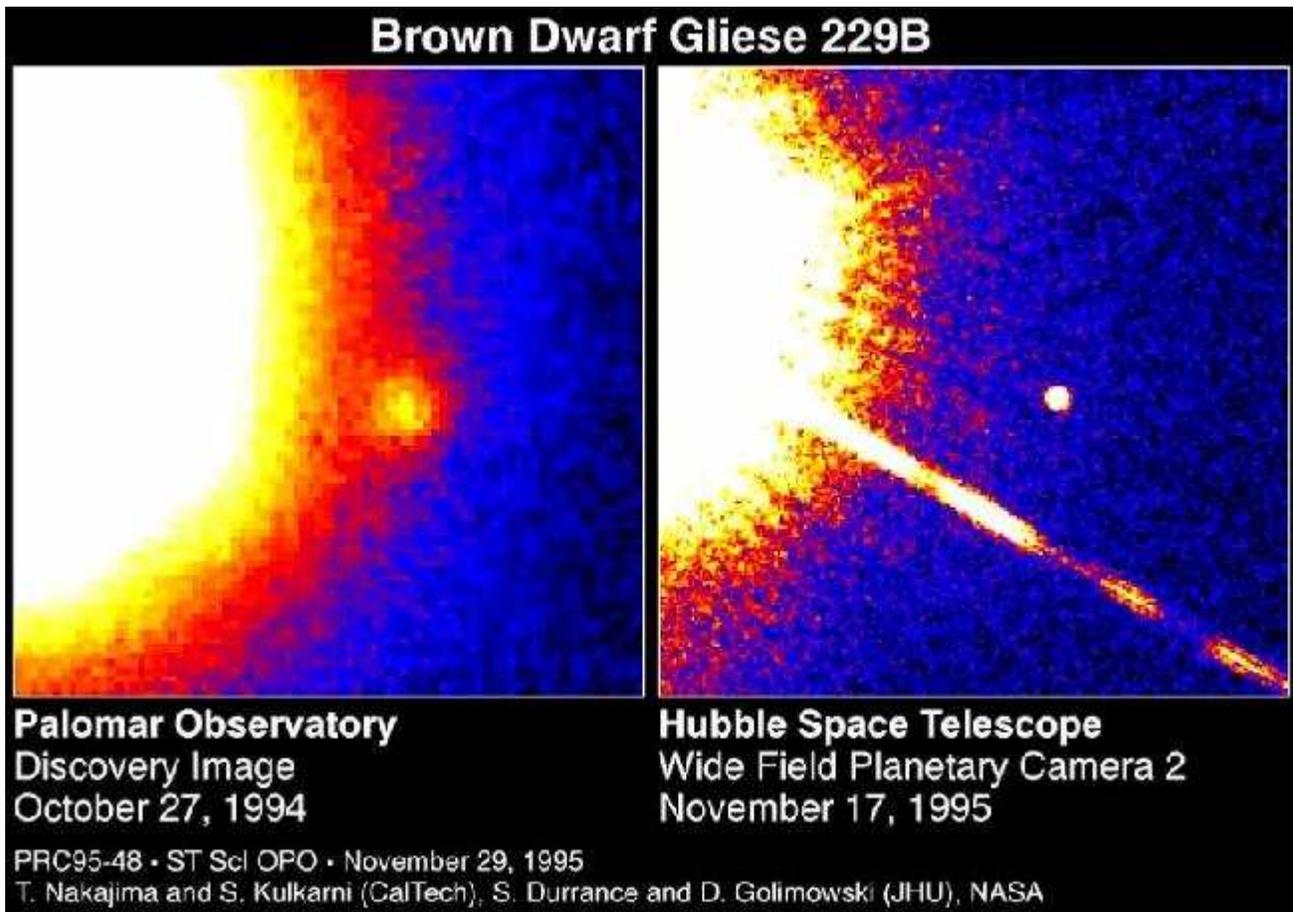
HST imaging found that Betelgeuse is one of the largest stars, almost the size of our whole solar system.



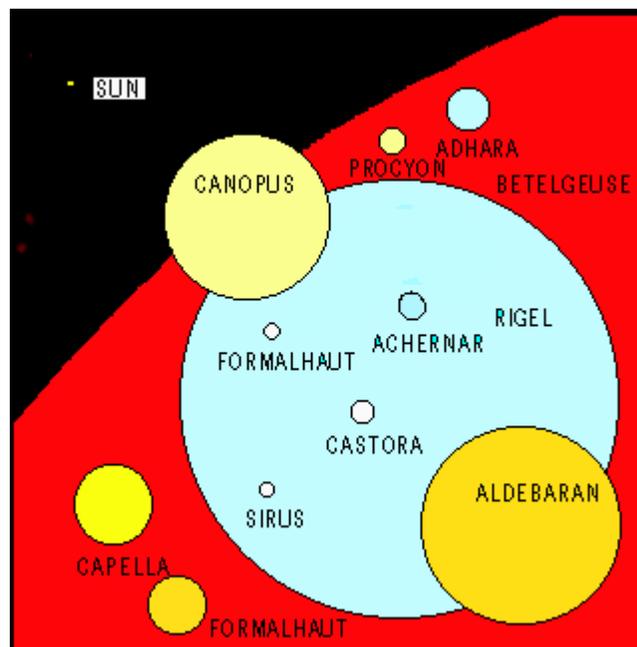
The other extreme was also found, that there exist a class of very small stars called white and brown dwarfs, with sizes close to the size of the Earth. White dwarfs are faint, but hot (=blue).



Brown dwarfs are very faint and quite red (=cool) making them hard to find in the sky.



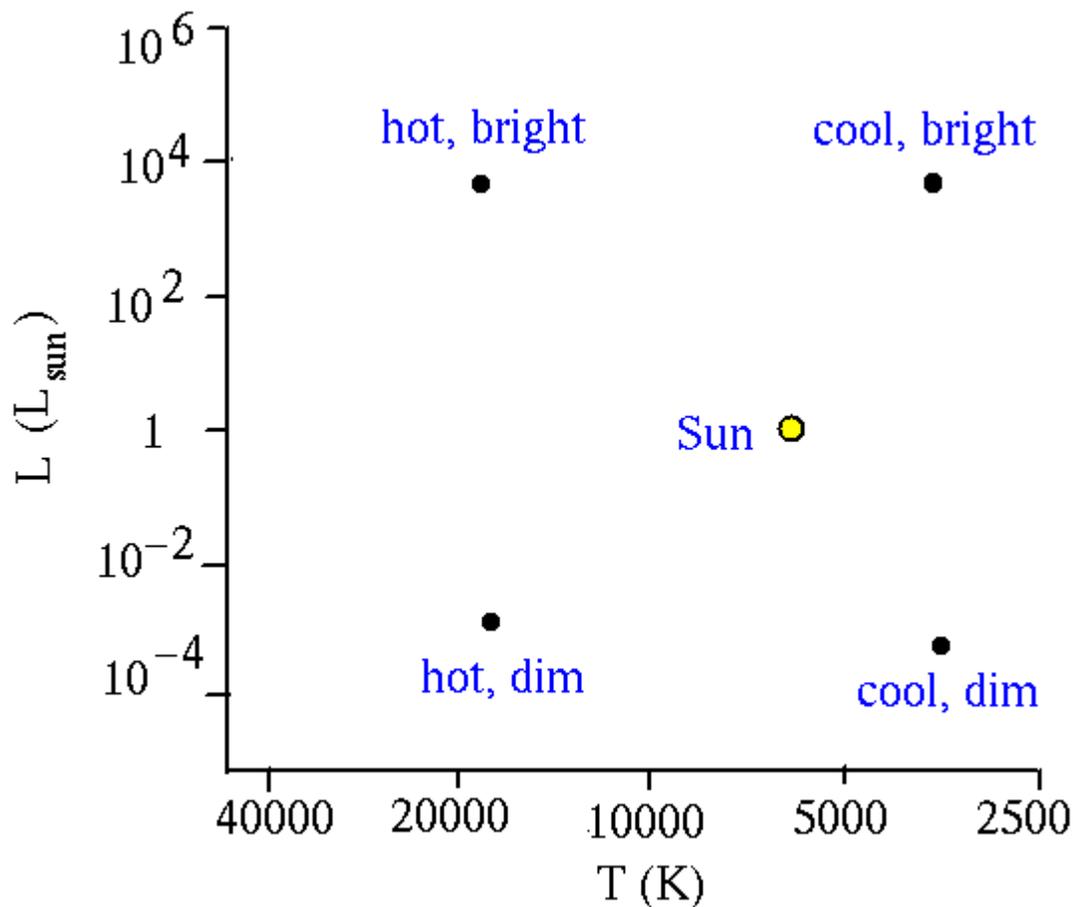
Red and blue supergiant stars, as well as giant stars exist. The following is a comparison of these types.



h) The Hertzsprung-Russell Diagram

In 1905, Danish astronomer Einar Hertzsprung, and independently American astronomer Henry Norris Russell, noticed that the luminosity of stars decreased from spectral type O to M. They developed the technique of plotting absolute magnitude for a star versus its spectral type to look for families of stellar type.

These diagrams, called the Hertzsprung-Russell or HR diagrams, plot luminosity in solar units on the Y axis and stellar temperature on the X axis, as shown below.

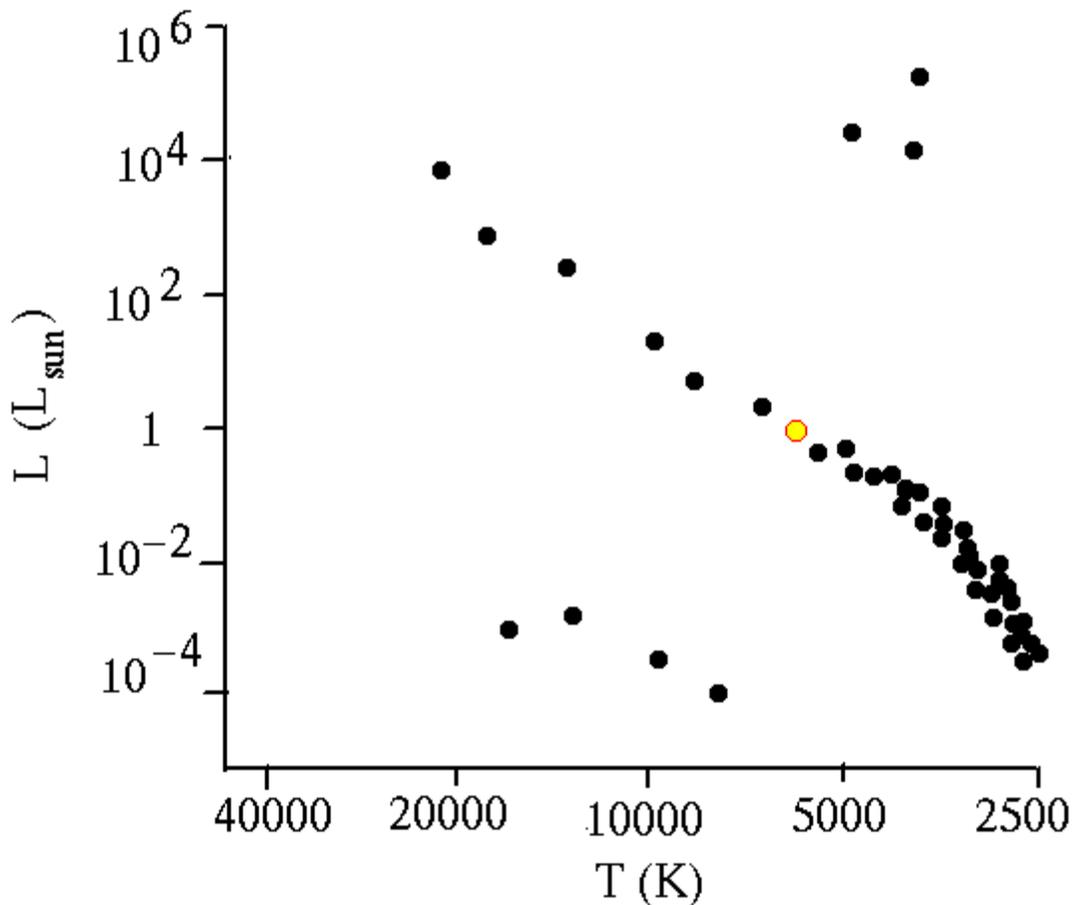


Notice that the scales are not linear. Hot stars inhabit the left hand side of the diagram, cool stars the right hand side. Bright stars at the top, faint stars at the bottom. Our Sun is a fairly average star and sits near the middle.

Notice that the vertical scale is luminosity as fractions of a stars absolute luminosity compared to the Sun. A star that is similar in brightness to the Sun has a L value of 1. A star that is 10,000 times brighter than the Sun has a L value of 10⁴. One that is 100 times fainter than the Sun has a L value of 10⁻². Likewise the horizontal axis is plotted in terms of temperature as measured in Kelvins. By custom we place hot stars on the left

and cool stars on the right, backwards from a normal graph.

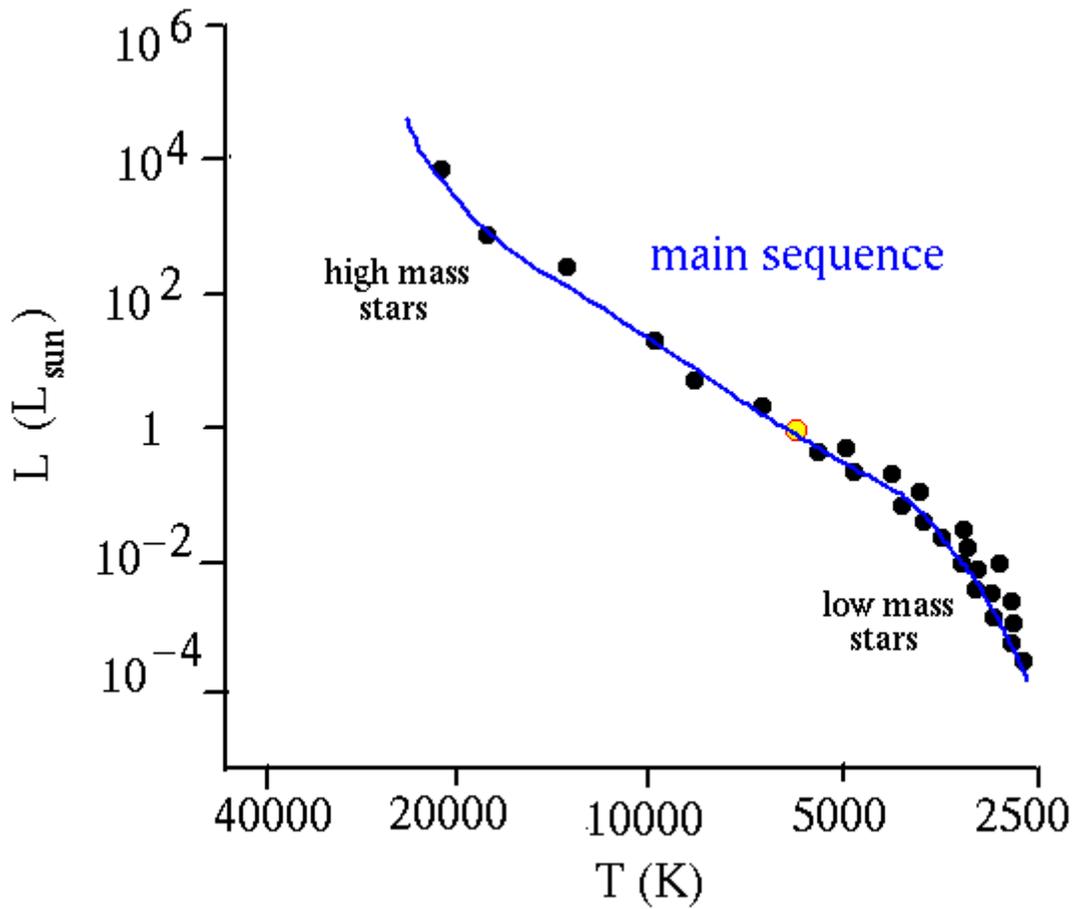
The HR diagram became one of our most powerful tools for understanding the physics of stars. Note, however, that an astronomer needs to measure two colors from the light of a particular star to get its temperature, then you need to know the distance to the star by parallax to convert its apparent magnitude into its absolute magnitude. Thus, the study of stars was a complicated and detailed process in the early days. The simplest stars to understand were the ones near our Solar System, with the largest parallaxes and the brightest apparent magnitudes. A plot of the nearest stars on the HR diagram is shown below:



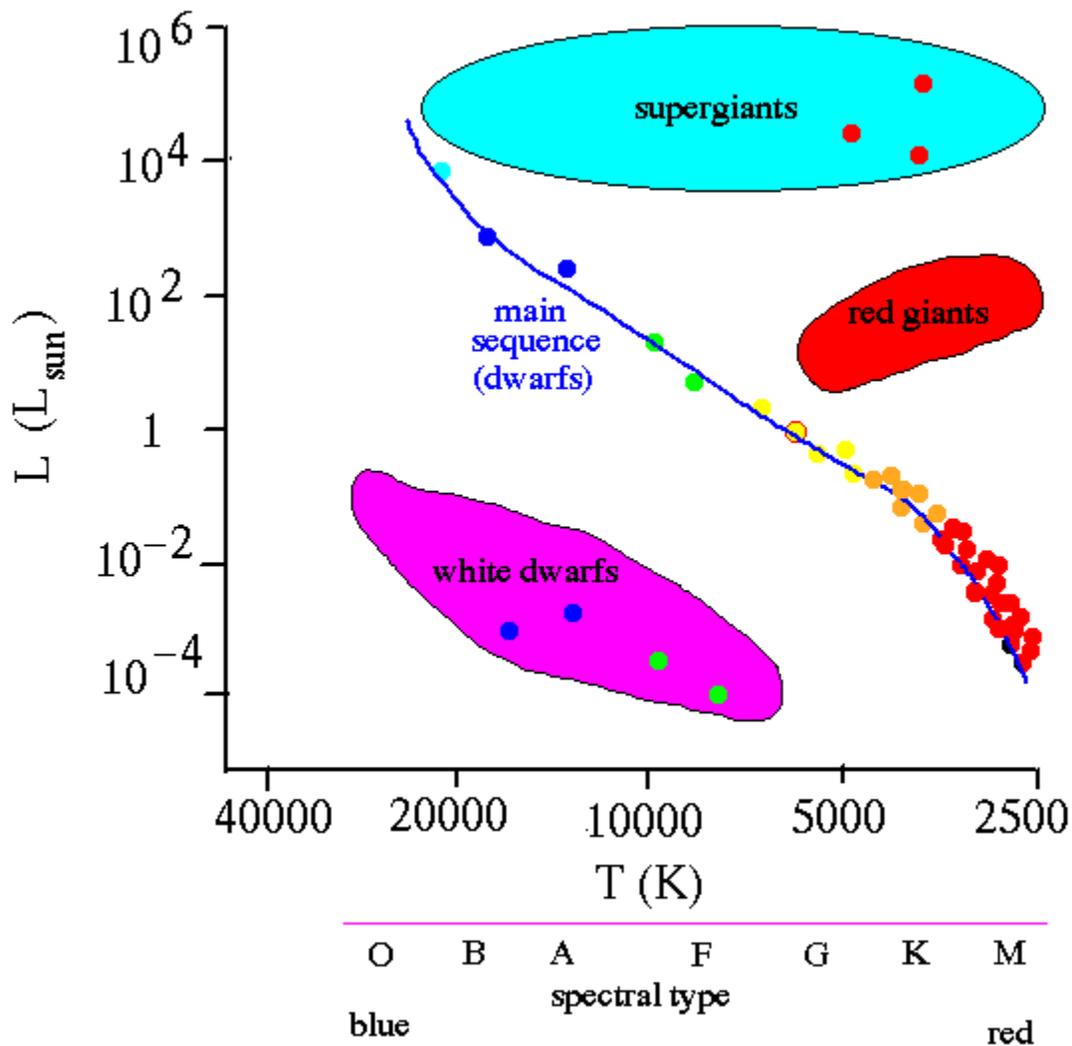
Most stars in the solar neighborhood are fainter and cooler than the Sun. There are also a handful of stars which are red and very bright (called red supergiants) and a few stars that are hot, but very faint (called white dwarfs). We will see in a later lecture that stars begin their life on the main sequence then evolve to different parts of the HR diagram.

Most of the stars in the above diagram fall on a curve that we call the main sequence. This is a region where most normal stars occur. Normal, in astronomy terms, means that

they are young (a few billion years old) and burning hydrogen in their cores. As time goes on, star change or evolve as the physics in their cores change. But for most of the lifetime of a star it sits somewhere on the main sequence.



Several regions of the HR diagram have been given names, although stars can occupy any portion. The brightest stars are called supergiants. Star clusters are rich in stars just off the main sequence called red giants. Main sequence stars are called dwarfs. And the faint, hot stars are called white dwarfs.

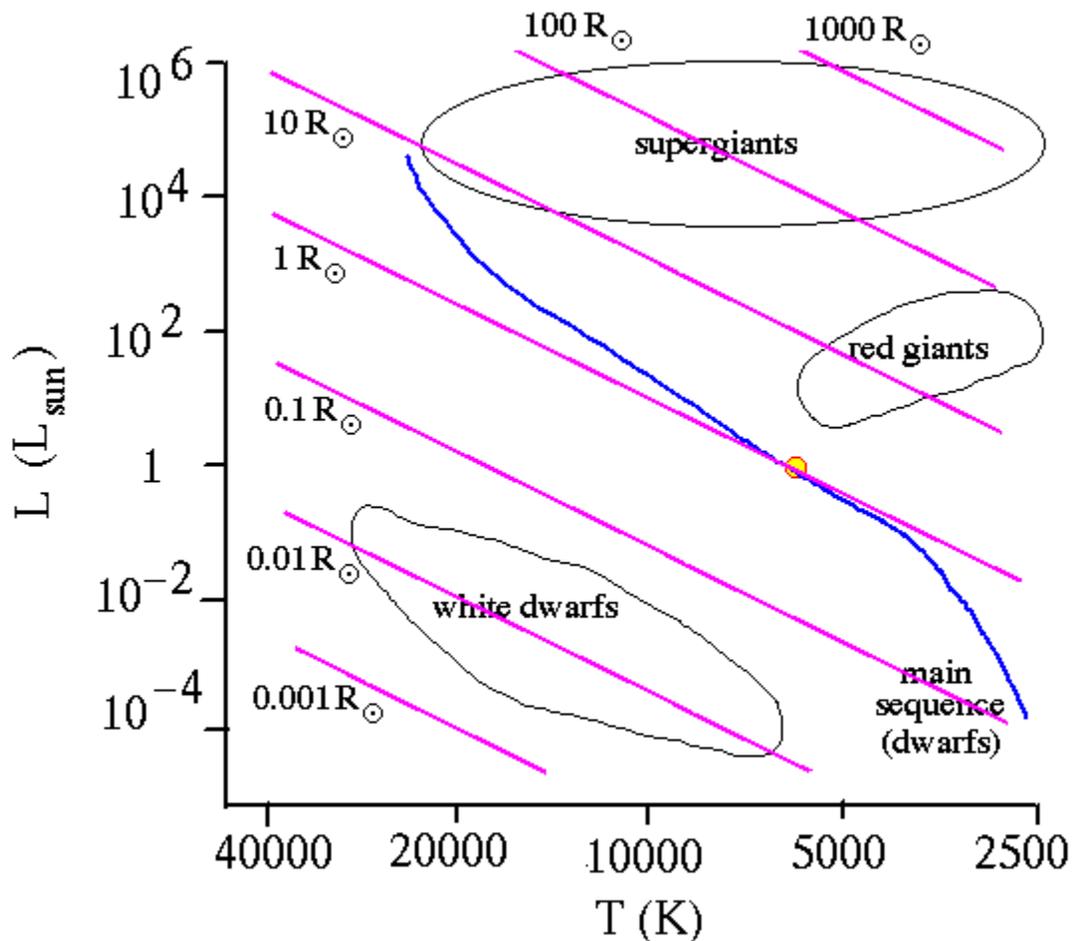


The spectral classification types were more accurate than attempts to measure the temperature of a star by its color. So often the temperature scale on the horizontal axis is replaced by spectral types, OBAFGKM. This had the advantage of being more linear than temperature (nicely spaced letters) and contained more information about the star than just its temperature (the state of its atoms).

The HR diagram becomes a calculating tool when one realizes that temperature, luminosity and size (radius) are all related by Stefan-Boltzmann's law.

Knowing from laboratory measurements that Stefan-Boltzmann's constant is 5.67×10^{-8} allows one to calculate the luminosity of a star in units of watts (like a light bulb) if we know the radius of the star in meters and the temperature in kelvins. For example, the Sun is 6.96×10^8 meters in radius and has a surface temperature of 5780K. Therefore, its luminosity is 3.84×10^{26} watts.

On a log-log plot, the R^2 term in the above equations is a straight line on an HR diagram. This means that on a HR diagram, a star's size is easy to read off once its luminosity and color are known.

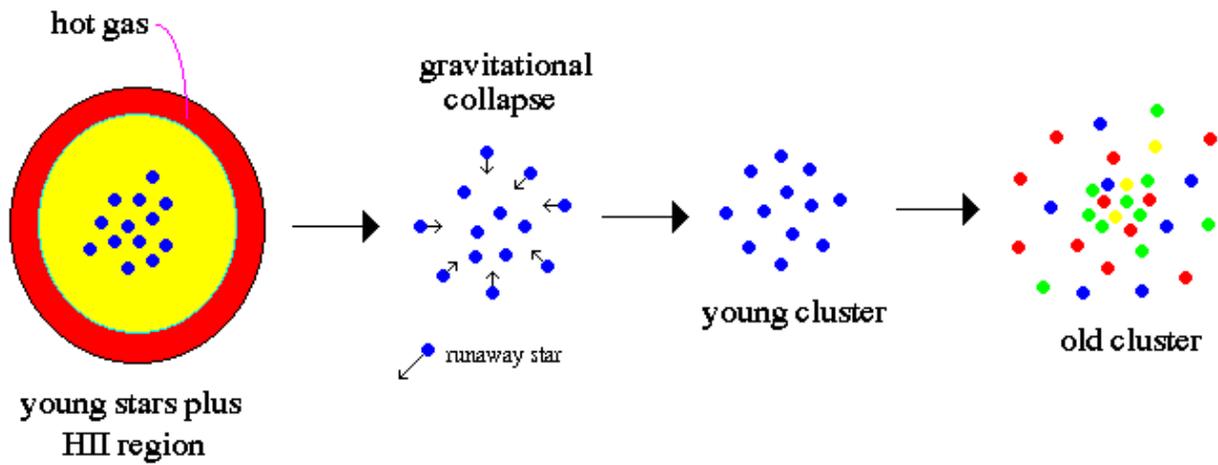


The HR diagram is a key tool in tracing the evolution of stars. Stars begin their life on the main sequence, but then evolve off into red giant phase and supergiant phase before dying as white dwarfs or some more violent endpoint.

i) Star Clusters

When stars are born they develop from large clouds of molecular gas. This means that they form in groups or clusters, since molecular clouds are composed of hundreds of solar masses of material. After the remnant gas is heated and blown away, the stars collect together by gravity. During the exchange of energy between the stars, some stars reach escape velocity from the protocluster and become runaway stars. The rest become

gravitationally bound, meaning they will exist as collection orbiting each other forever.



When a cluster is young, the brightest members are O, B and A stars. Young clusters in our Galaxy are called open clusters due to their loose appearance. They usually contain between 100 and 1,000 members. One example is the binary cluster below:



And the Jewel Box cluster:



Early in the formation of our Galaxy, very large, globular clusters formed from giant molecular clouds. Each contain over 10,000 members, appear very compact and have the oldest stars in the Universe. One example is M13 (the 13th object in the Messier catalogs) shown below:

Globular Cluster NGC 6093



Hubble
Heritage

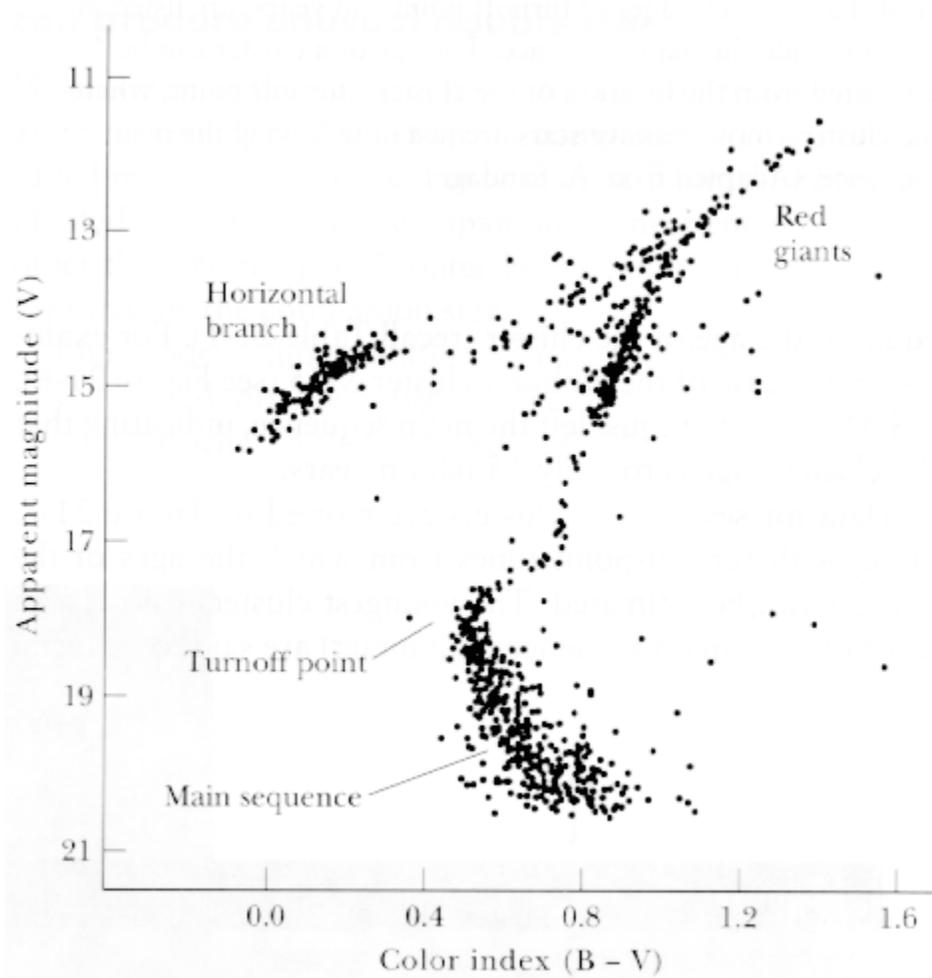
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Since all the stars in a cluster formed at the same time, they are all the same age. A very young cluster will have a HR diagram with a cluster of T-Tauri stars evolving towards the main sequence. As time passes the most massive stars at the top of the main sequence

evolve into red giants. Therefore, the older the cluster, the fewer stars to be found at the top of the main sequence, and an obvious grouping of red giants will be seen at the top right of the HR diagram.

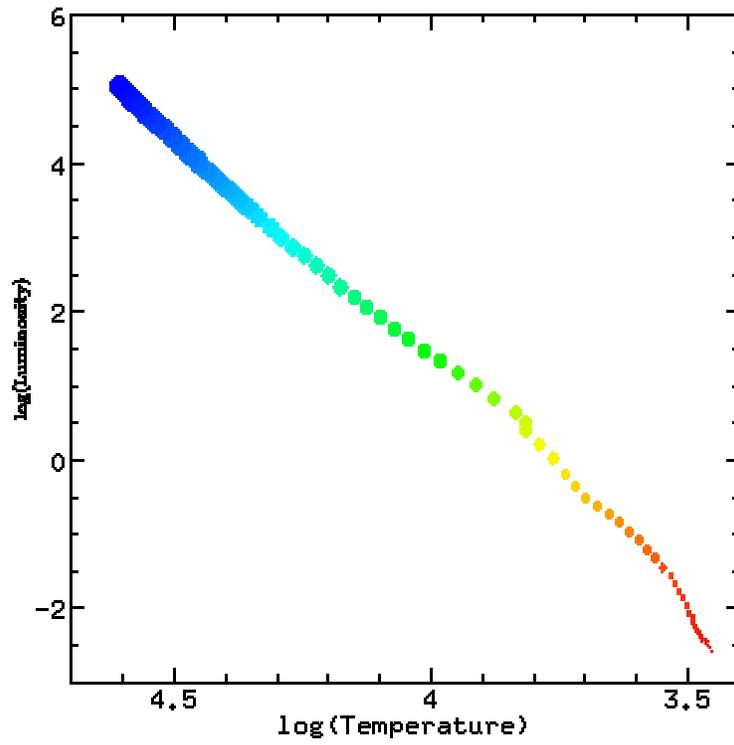
This effect, of an evolving HR diagram with age, becomes a powerful test of our stellar evolution models. A computer model can be built that follows the changes in stars of various masses with time. Then an theoretical HR diagram can be built at each timestep. Observations of star clusters are compared to these computer generated HR diagrams to test of our understanding of stellar physics.

Observations of star clusters consist of performing photometry on as many individual stars that can be measured in a cluster. Each star is plotted by its color and magnitude on the HR diagram. Shown below is one such diagram for the globular cluster M13.

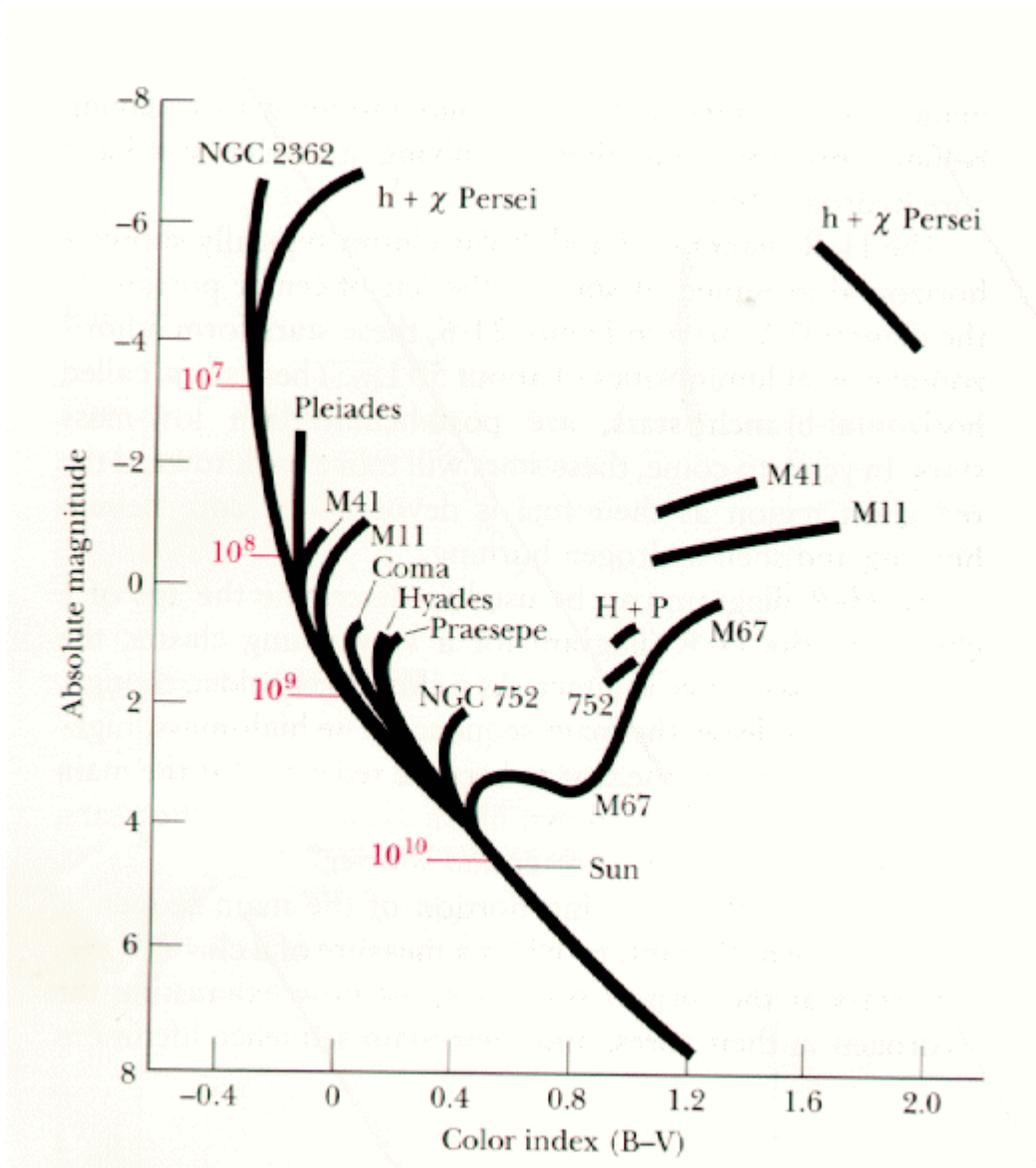


Note that the main sequence only exists for low mass G, K and M stars. All stars bluer than the turn-off point have exhausted their hydrogen fuel and evolved into red giants millions and billions of years ago. Also visible is a clear red giant branch and a post-red

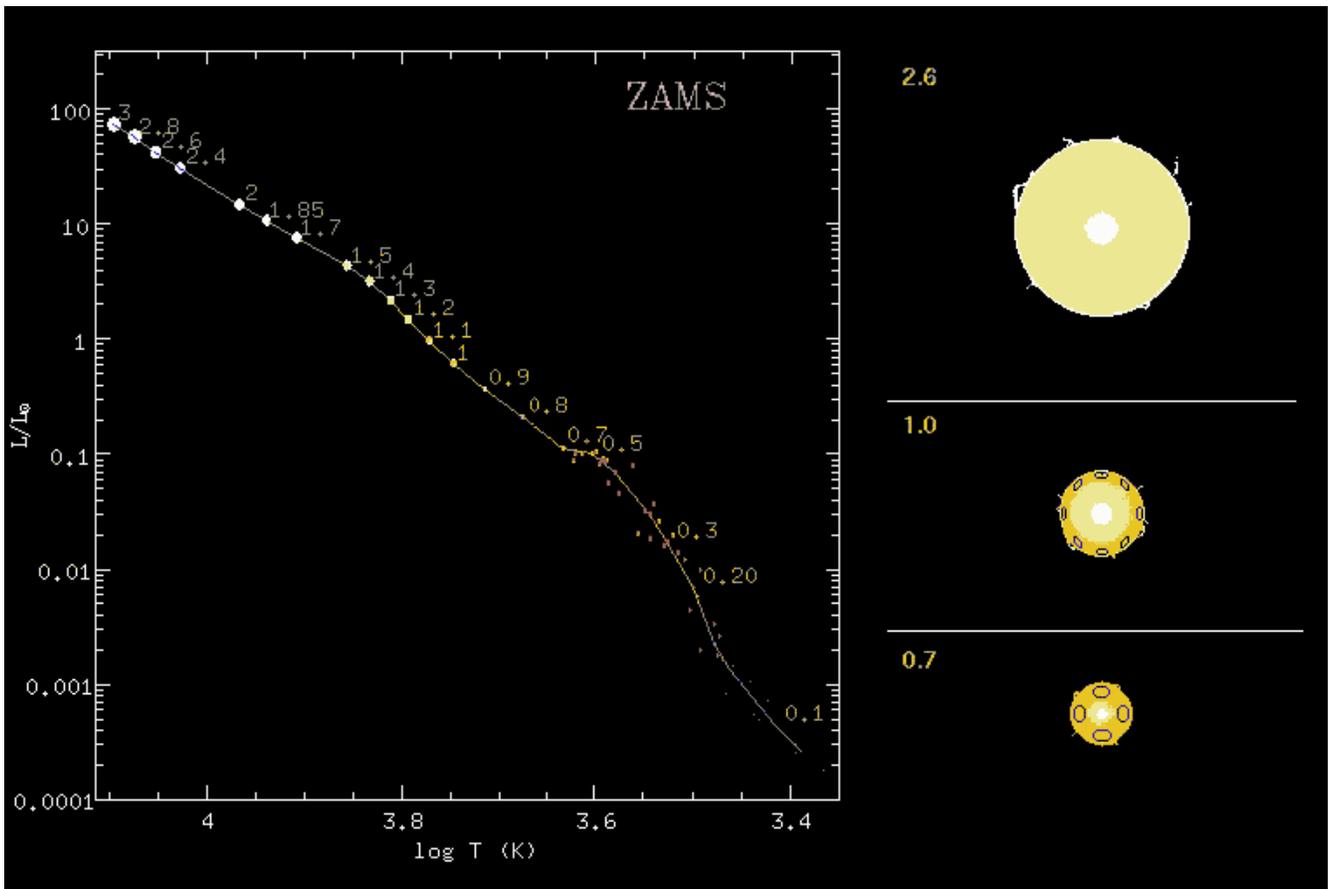
giant phase region, the horizontal branch.



Plotting various star cluster HR diagrams together gives the following plot



Understanding the changes in the lifetime of a main sequence star is a simple matter of nuclear physics, where we can calibrate the turn-off points for various clusters to give their ages. This, then, provides a tool to understand how our Galaxy formed, by mapping the positions and characteristics of star clusters with known ages. When this is done it is found that old clusters form a halo around our Galaxy, young clusters are found in the arms of our spiral galaxy near regions of gas and dust.



The above animation is a detailed computer simulated HR diagram. The horizontal axis is the log of temperature, the y axis is the luminosity of the stars. The main sequence is the line through the middle of the figure. Notice the stars evolving off the main sequence labeled by their masses in solar masses, the high mass ones first. To the right is a diagram of the interiors of three stars, masses 2.6, 1.0 and 0.7 solar masses. The cutaways shows what is happening inside the stars as they evolve. Notice the enlarging envelopes as the stars become red giants.